Non-equilibrium helium ionization in the solar atmosphere

by

Thomas Peter Golding

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Institute of Theoretical Astrophysics
Faculty of Mathematics and Natural Science
University of Oslo
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Part I

Background
The solar atmosphere is, roughly speaking, the part of the sun from which emitted light is able to escape. It is traditionally split into four different regions: the photosphere, the chromosphere, the transition region, and the corona. Figure 1.1 shows the temperature and the mass density of a column taken from a numerical model of the solar atmosphere (see Section 2.2 for information about the model). The vertical extent of the four atmospheric regions is indicated in the figure. The figure shows the conditions from the chosen atmospheric column at a particular time. At a later time or in a

![Figure 1.1: Model of the solar atmosphere showing the temperature (solid line, scales to the left) and the mass density (dashed line, scales to the right). The data is taken from a column of a numerical simulation.](image)
Figure 1.2: The solar atmosphere shown in two spectral positions in the H-α line. The left panel shows the atmosphere in the far blue wing at -60 km/s from line center which is shown in the right panel. These observations were made on the 15th of June 2008 by the CRISP instrument at the SST.

different column, the extent of the four atmospheric regions will be different. The terms photosphere, chromosphere, transition region, and corona are therefore acting more as indicators of particular physical regimes rather than indicators of particular heights in the atmosphere. The next sections contain a brief description of the different regimes of the atmosphere.

The photosphere

The solar photosphere is the lower part of the solar atmosphere. It is roughly 500 km thick. In the layers below, energy is transported upwards by convective motions in the convection zone. The photosphere is where the convection zone gas emerges, cools off radiatively, and sinks back down. The visible pattern of the convection process is known as granulation. The left panel of Figure 1.2 shows the granular pattern in an observation of the H-α line done with the CRISP instrument (Scharmer et al., 2008) mounted on the Swedish 1-m Solar Telescope (SST) (Scharmer et al., 2003). The bright points in the granulation correspond to locations with strong magnetic field concentrations sticking out of the photosphere (Dunn & Zirker, 1973; Stein, 2012). Outside of these bright patches the magnetic field strength is lower. In the photosphere gas pressure forces are for the most part larger than the magnetic forces. This means that the gas pressure dominates the dynamics of the photosphere. The temperature of the photosphere is around 5000 K.
At such temperatures the gas is mostly neutral. The radiative cooling of the photospheric gas happens because the opacity of the material at the visual light wavelengths becomes low enough for photons to escape. This marks the transition where the radiation temperature decouples from the gas temperature. Below this point, where the density is high and the radiation temperature equals the gas temperature, the ionization and excitation state can be described by local thermodynamic equilibrium (LTE).

The chromosphere

The chromosphere is the layer above the photosphere. Here the gas has a lower density by orders of magnitude, resulting in a lower gas pressure. The relative importance of the magnetic forces therefore increases and the magnetic forces can dominate the dynamics. An example of a magnetic feature are chromospheric fibrils, shown in the right panel of Figure 1.2, who are correlated with the magnetic field (de la Cruz Rodríguez & Socas-Navarro, 2011; Leenaarts et al., 2015).

The chromosphere is a dynamic region. Waves are excited in deeper layers and propagate upwards (Carlsson & Stein, 1997; Wedemeyer et al., 2004; Morton et al., 2012). Chromospheric jets known as spicules (Beckers, 1968) are occurring everywhere all the time. Such dynamic features affect the thermodynamic state of the chromosphere. The thermodynamic state is closely related to the ionization state of the gas. The low density results in low collisional transition rates, and the radiative transitions become important. The lower transition rates lead to ionization-recombination timescales comparable to the dynamical timescales. Therefore the ionization state of the chromosphere is not in equilibrium (Carlsson & Stein, 2002; Leenaarts et al., 2007).

The transition region

The transition region is a thin layer of the atmosphere where the temperature increases from around $10^4$ K to $10^6$ K. Moving upward from the chromosphere the gas is heated, and much of this energy goes in to ionizing the atoms. When all atoms are ionized the gas can no longer deposit energy in ions and the temperature rises. Figure 1.3 shows an Interface Region Imaging Spectrograph (IRIS) 1400 slit-jaw image (De Pontieu et al., 2014). The image gives an impression of what the transition region looks like since it is dominated by lines forming at around 60,000 kelvin.

The corona

The corona is the upper part of the atmosphere. Here the temperature is of the order of $10^6$ K. It is highly ionized, and the magnetic forces completely dominate the dynamics. How the chromosphere and corona is heated to
Figure 1.3: Slit-jaw image 1400 from IRIS, taken on the 6th of November 2014. The image shows the transition region. Si IV lines forming at a temperature roughly 60,000 K are dominating in the 1400 passband.

maintain such high temperatures is a topic of debate, but magnetic fields are believed to be central (Klimchuk, 2006).
Modelling the solar atmosphere

Information about the solar atmosphere is carried by the light it emits. Imprinted in the spectrum lie clues about the physical conditions under which the light was formed. The formation of the solar spectrum is complex. Therefore, using measurements of a solar spectrum to diagnose the physical conditions is a challenge. For this purpose, models of the solar atmosphere are of great value. The work presented in this thesis was carried out using different computer codes to model properties of the solar atmosphere. This chapter contains the basic equations used in these codes, and it points out problems related to their solution.

2.1 Physical model

The radiation-magnetohydrodynamics equations

Following a fluid approach, the equations governing the solar atmosphere are the equations of radiation-magnetohydrodynamics (RMHD). They are expressed,

\[ \frac{\partial \rho}{\partial t} = - \nabla \cdot (\rho u) \]  
\[ \frac{\partial (\rho u)}{\partial t} = - \nabla \cdot (\rho uu - \epsilon) - \nabla P + J \times B + \rho g \]  
\[ \mu J = \nabla \times B \]  
\[ \frac{\partial B}{\partial t} = \nabla \times (u \times B - \eta J) \]  
\[ \frac{\partial e}{\partial t} = - \nabla \cdot (eu) - P \nabla \cdot u + Q_{\text{rad}} + Q_{\text{other}} \]

where \( \rho, u, \epsilon, P, J, B, g, \mu, \eta, e, \) and \( Q_{\text{rad}} \) are the mass density, flow velocity, stress tensor, gas pressure, current density, magnetic field strength, magnetic field intensity, and energy generation, respectively.
gravitational acceleration, vacuum permeability, magnetic diffusivity, internal energy, and radiative heating rate, respectively. $Q_{\text{other}}$ is the heating rate due to heat conduction, viscous dissipation, and ohmic dissipation.

### Equation of state

Additional equations relate the mass density and internal energy. These are

\[
P = \left( \sum_{i,k} n_{ik} + n_e \right) k_B T \tag{2.6}
\]

\[
e = \frac{3}{2} \left( \sum_{i,k} n_{ik} + n_e \right) k_B T + \sum_{i,j} n_{ik} E_{ik} \tag{2.7}
\]

\[
n_e = \sum_{ik} n_{ik} Y_{ik} \tag{2.8}
\]

\[
\sum_i n_{ik} = A_k \rho. \tag{2.9}
\]

Here the index $k$ indicates the element and $i$ indicates the atomic state, so $n_{ik}$ and $E_{ik}$ denote the population density and energy of an atom of element $k$ in state $i$. $Y_{ik}$ is the number of electrons lost from an atom of type $k$ in state $i$ compared to a neutral state. $n_e$ is the electron density, $T$ is the temperature, and $k_B$ is the Boltzmann constant.

The pressure, temperature and electron density can be computed once the population densities, $n_{ik}$, are determined. In general these densities are solutions to a set of rate equations:

\[
\frac{\partial n_{ik}}{\partial t} = -\nabla \cdot (u n_{ik}) + \sum_{j \neq i} n_{jk} P_{ijk} - n_{ik} \sum_{j \neq i} P_{ijk}. \tag{2.10}
\]

Here $P_{ijk}$ is the probability of an atom of type $k$ in the state $i$ transitioning to the state $j$, given by the sum of a collisional ($C_{ijk}$) and radiative ($R_{ijk}$) transition coefficient. These are differential equations that should be integrated forward in time together with the RMHD equations, making the population densities at any given time dependent on the history of the atmosphere.

### Radiative transfer

Radiative transfer is needed to obtain the radiative heating and solve the rate equations. The radiative heating is given by the frequency integral

\[
Q_{\text{rad}} = 4\pi \int_0^\infty \chi_\nu (J_\nu - S_\nu) \, d\nu, \tag{2.11}
\]
where $\chi_\nu$, $J_\nu$, and $S_\nu$ are the frequency dependent opacity, mean intensity, and source function. The radiative transition coefficient is given by a frequency integral involving the mean intensity,

$$R_{ijk} = \int_0^\infty J_\nu f(\nu) \, d\nu,$$

(2.12)

where the function $f(\nu)$ is dependent on the which atom is under consideration and different for bound-bound and bound-free transitions. $J_\nu$ is the angular average of the specific intensity, $I_\nu$, which is the solution of the equation of radiative transfer,

$$\frac{dI_\nu}{d\tau_\nu} = S_\nu - I_\nu,$$

(2.13)

where $\tau_\nu$ is the frequency dependent optical path length along a ray.

### 2.2 Bifrost, a numerical model

The stellar atmosphere code Bifrost (Gudiksen et al., 2011) is central to the work presented in this thesis, and a brief description of the code is given here. The code solves the RMHD equations (2.1-2.5) in parallel on a staggered cartesian grid in two or three spatial dimensions. The computational domain is split up in sub-domains which are distributed among the running CPUs. The heat conduction follows Spitzer (1956) and is oriented along the magnetic field lines. The radiative transport needed for the calculation of the radiative heating rate, $Q_{rad}$, is done in three different modules of the code. The first, dealing with the radiation transport in the convection zone, photosphere and low chromosphere, is utilizing an opacity binning method (Nordlund, 1982), including scattering (Skartlien, 2000; Hayek et al., 2010). The second, dealing with losses from strong chromospheric lines, is approximated by using various fitting functions derived from 1D hydrodynamical simulations (Carlsson & Leenaarts, 2012). The third, dealing with the cooling from EUV lines in the thin transition region and corona, is found assuming coronal ionization equilibrium, in which case the cooling function can be determined by the software package CHIANTI (Dere et al., 1997; Del Zanna et al., 2015).

### Equation of state

The pressure, temperature and electron density are determined by equations 2.6-2.9, which depend on the atomic number densities. The general approach is to determine the atomic number densities by solving a set of rate equations (2.10). This requires knowledge of the transition rate coefficients. While the contributions due to collisional transitions are expressed as functions of
temperature and electron density (in addition to atomic data), the radiative transition coefficients pose a problem as they are given by frequency integrals over the mean intensity weighted by some function (eq. 2.12).

The rate equations (2.10) express that any change in atomic number densities are caused by either advection (\( \nabla \cdot \mathbf{u} n_{ik} \)) or atoms making transitions from one state, \( i \), to a different state, \( j \), (\( n_{ik} P_{ijk} \)). In the convection zone and photosphere the rate of change due to atomic transitions is faster than the rate of change due to advection. The gas is optically thick, which couples the radiation to the local temperature. Under such conditions the collisional and radiative rate coefficients ensure instant equilibration to the LTE ionization and excitation state. In this case the atomic number densities can be computed from the simpler Saha-Boltzmann equations (Mihalas, 1978). In the chromosphere and above, the gas becomes increasingly transparent for photons and non-local coupling through radiation becomes important. Also, the transition rate coefficients decrease in value. This increases the timescale of atomic transitions. Therefore LTE is not valid, and the full solution of the set of non-equilibrium rate equations is required.

Modern numerical codes, such as Bifrost, running on contemporary computer facilities are not able to solve the RMHD equations including a detailed account of multidimensional multilevel radiative transfer, resolving individual line profiles. This has prevented inclusion of non-equilibrium ionization in such simulations, and assuming LTE has been the the common way of approximating the equation of state (e.g. Freytag et al., 2012; Hansteen et al., 2015; Wray et al., 2015; Martínez-Sykora et al., 2016). However, the calculation of non-equilibrium atomic number densities needed for a general solution of the RMHD equation set can be simplified without a too large loss of realism. These simplifications are an important theme of this thesis.

**Non-equilibrium ionization**

Taking non-equilibrium ionization of abundant elements into account when modelling the solar atmosphere has a significant impact on the energy balance of the gas. The most abundant elements in the solar gas composition are hydrogen and helium (Grevesse & Sauval, 1998), and they dominate the equation of state. Consider neutral solar gas at temperature 10000 K. It consists of 90% hydrogen and 10% helium. Doubling the temperature, ionizing 7% of the hydrogen, or ionizing 35% of the helium, all require the same amount of energy. The ionization state of other less abundant elements have a negligible effect on the energy balance.

**Hydrogen**

Non-equilibrium ionization of hydrogen results in hotter shock fronts and colder post shock gas in the chromosphere due to long equilibration timescales
2.3 Calculating the emergent intensity

(Carlsson & Stein, 1992, 2002). Bifrost has an option for including non-equilibrium ionization of hydrogen (Leenaarts et al., 2007, 2011). When the option is switched on, the code solves the rate equations for five neutral states, the ionized state, and one molecular state of hydrogen. The rate equations are solved together with conservation equations for internal energy (2.7), charge (2.8), and hydrogen particles (2.9). The method exploits prescribed radiative transition rate coefficients (Sollum, 1999).

Helium

Helium is mostly neutral in the photosphere, partially ionized in the chromosphere and transition region, and fully ionized in the corona. Describing the ionization state of helium and the effects on the solar atmosphere is a central theme of this PhD project and is elaborated in chapter 4.2.

2.3 Calculating the emergent intensity

Measurements of the intensity emerging from the solar atmosphere are the primary quantities for studies of the solar atmosphere. As already mentioned, Bifrost only computes the intensity for a few frequency bins. To be able to directly compare individual line profiles and continua, a resolution of a few hundred frequency points are needed. The procedure of computing the emergent intensity is done by a separate compute code. In this work the code Multi3d (Leenaarts & Carlsson, 2009) is used. Multi3d can solve the non-LTE radiative transfer problem on a single snapshot from a Bifrost simulation. That means it will solve the radiative transfer equation together with the rate equations. Statistical equilibrium (SE) is assumed, which means that the terms $\frac{\partial n_{ik}}{\partial t}$ and $\nabla \cdot (u n_{ik})$ are both set to zero.

There are, however, situations where the SE assumption is not fully justified. Possible non-equilibrium effects will be missed in situations where the desired diagnostic is formed in a region where the rates are slow and the dynamics are fast. A method of approximating non-equilibrium non-LTE radiative transfer was developed as part of this thesis and a brief description is given in section 4.3.
Helium diagnostics

Helium spectral lines are popular diagnostics of the solar atmosphere. Helium resonance EUV lines, which are not observable from the ground, have been, or are currently, observed by SOHO/CDS (Harrison et al., 1995), SOHO/SUMER (Wilhelm et al., 1995), HINODE/EIS (Culhane et al., 2007), and SDO/AIA (Lemen et al., 2012). The number one helium line observed from the ground is the $\lambda 10830$ of He I. Examples of instruments in operation observing this line are VTT/TIP II (Collados et al., 2007), NST/NIRIS (Cao et al., 2012), and GREGOR/GRIS (Collados et al., 2012).

3.1 Helium EUV lines

The formation of the helium EUV lines have been debated for some decades. Jordan (1975) pointed out that the observed line intensities of the first resonance lines of He I and He II were higher by a factor of $\sim 10$, compared to what her emission measure model predicted. The same trend has been confirmed in other studies (MacPherson & Jordan, 1999; Pietarila & Judge, 2004; Giunta et al., 2015). Broadly speaking there are two mechanisms that have been proposed to explain the high EUV line intensities. The first one was proposed already by Jordan (1975). She suggested a mechanism admitting cold atoms and ions to be excited by hot electrons. This would enhance the intensities of the lines. Drivers of such a mechanism have been suggested to be microturbulent motions Jordan (1980); Andretta et al. (2000); Smith & Jordan (2002); Pietarila & Judge (2004), particle diffusion Fontenla et al. (1990, 1991, 1993), fluid flows Fontenla et al. (2002), and non-maxwellian electron distributions Smith (2003). All of these mechanism drivers are able to enhance the intensity of the helium EUV lines. The most popular one seems to be the one related to microturbulent fluid motions. This is often referred to as velocity redistribution or VR for short, after Andretta et al. (2000). The other mechanism that has been proposed is
the so called photoionization-recombination or PR mechanism (Zirin, 1975, 1988, 1996). Under a PR scenario a helium atom is photoionized, after which it recombines and cascades down to the ground state. The final step in the cascade will produce an EUV photon. Theoretical investigation of this mechanism have shown it to enhance the He I \( \lambda 584 \) line under quiet conditions, but less so in more active regions (Andretta & Jones, 1997; Mauas et al., 2005).

### 3.2 He I \( \lambda 10830 \)

The He I \( \lambda 10830 \) spectral line is perhaps the most commonly observed helium line from ground-based observatories. Theoretical modelling efforts have shown that the line forms in the chromosphere and that it is sensitive to the incident radiation from the transition region and corona (Fontenla et al., 1993; Avrett et al., 1994; Andretta & Jones, 1997; Mauas et al., 2005; Centeno et al., 2008). In other words it carries information not only about chromospheric conditions, but also about the activity of the regions where the EUV photons are emitted. In addition, the line is sensitive to both the Hanle and Zeeman effect, and is as such a good tool for measuring the chromospheric magnetic field (Rueedi et al., 1995; Wiegelmann et al., 2014; Lagg et al., 2015).

### 3.3 Modern models are needed

The modelling of helium lines emerging from the solar atmosphere has so far been based on assumptions of static or stationary conditions as well as simple geometry. Observational facilities are getting better and better. While comparing modern observations to 1D static models is good, supplementing with 3D time dependent models is better. A part of this PhD project is to create such modern models, and this is elaborated in section 4.3.
CHAPTER 4

PhD project summary

4.1 The project goals

The goal of the PhD project presented in this thesis was to describe how non-equilibrium helium ionization affects the energetics and helium line formation in the context of the solar atmosphere. First a computer code was developed, capable of producing numerical models of the solar atmosphere where non-equilibrium ionization of abundant elements was taken into account. From such models the effects on the energy state was studied. These models also served as a suitable basis for studies of helium line formation with the help of radiative transfer codes. The rest of this chapter is devoted to the work that was done. Section 4.2 gives a description of how helium ionizes in the solar atmosphere, what the effects are, as well as how this was implemented in Bifrost. Section 4.3 gives a summary of our findings regarding the formation of helium lines.

4.2 Helium ionization in solar atmosphere models

Helium ionization in a 1D model

The ionization state of helium is important for the energy balance of the solar atmosphere. Helium goes from mostly neutral in the photosphere to fully ionized in the corona. In the chromosphere and transition region helium is partially ionized. To assess which processes dominate in setting the ionization state of helium, we set up a solar atmosphere simulation using the 1D radiation-hydrodynamics code Radyn (Carlsson & Stein, 1992, 1995, 1997, 2002). The code simulates the physical model described in section 2.1, but without magnetic fields. Since Radyn solves in detail
the non-equilibrium rate equations (eq. 2.10), it is suitable for studies of ionization-recombination mechanisms. The simulation features waves excited in the lower boundary, steepening into shocks in the chromosphere. A constant incident EUV radiation field is imposed as the upper boundary condition for the specific intensity. The incident radiation represents the emission in EUV lines of the transition region and corona, and it is based on the models from Tobiska (1991). We use a 33 level helium model atom.

We find that the ionization state of helium in the chromosphere is set by photoionization due to the incident EUV radiation. In the transition region the temperature is high enough for collisions to contribute. The ionization is primarily counteracted by spontaneous radiative recombination to excited states, followed by a cascade process where the bound electron transitions down in a stepwise fashion ultimately reaching the ground state energy level \( n = 1 \). In the model the timescale of ionization-recombination in the chromosphere and transition region is comparable to the dynamic timescale of the atmosphere, leading to non-equilibrium helium ion fractions. The gas in the photosphere is dense and optically thick. The high density leads to high valued collisional transition rate coefficients. The optical thickness leads to a radiation field, and therefore also radiative transition rate coefficients, coupled to the local gas conditions. The high valued and locally determined transition rate coefficients result in LTE conditions.

**A simplified helium model atom**

A 3D RMHD code is unable to solve the multilevel non-LTE radiation transfer problem due to inadequate computing capacity available on contemporary systems. By reducing the complexity of the physical system, the problem can become tractable. We therefore construct a simplified 3 level helium model atom consisting of ground state He I, ground state He II, and He III. The role of all excited states is assumed to be temporary stepping stones in a recombination cascade. Based on the Radyn simulation we claim that this is a fair assumption. The idea is that all helium atoms that recombine to an excited state will de-excite down to the ground state, and they do not act as states from which collisional or radiative ionization is taking place. The role of the excited states can then be expressed by extra recombination rate from the He II and He III states. The recombination rate coefficients are tabulated and based on the values from the Radyn simulation. The 3 level model atom is simple enough to be used in a 3D RMHD code and complex enough to capture the essential physics. Details on the simplified model atom can be found in Golding et al. (2014) which is given in chapter 5.
The implementation of non-equilibrium helium ionization in Bifrost

We develop a method for including non-equilibrium helium ionization in Bifrost. The method exploits the simple 3 level helium model atom. The 3 atomic levels or states translate to 3 additional equations that must be solved. Two rate equations and one conservation equation for helium particles are chosen. These three equations are solved together with the nine equations of the non-equilibrium hydrogen package in the Bifrost code (see section 2.2). This way non-equilibrium hydrogen and helium ionization is taken into account simultaneously.

The EUV photons emitted from the transition region and corona, and their absorption in the helium and hydrogen continua, are self-consistently modelled. The emissivity is due to EUV lines from heavy ions. By assuming ionization equilibrium and optically thin conditions, given atomic data, the emissivity is expressed by a temperature dependent function multiplied by the hydrogen atom and electron densities. The opacity is taken to be a function of the number density of the ground states of He I, He II, and H I. The emissivity and opacity are computed for a few frequency bins only. Once the emissivity and opacity are known, the mean intensity field and transition rate coefficients can be calculated. Equipped with the radiation transfer quantities, the EUV part of the heating rate (eq. 2.11) and radiative rate coefficients for helium (eq. 2.12) are easily found.

Consequences of non-equilibrium ionization in Bifrost models

The non-equilibrium ionization equation of state removes a tendency of grid cells having certain preferred temperatures that are present when using an LTE equation of state. The grid cells in question are the ones corresponding to the chromosphere. The preferred temperature are manifested as horizontal dark bands in figure 4.1, which shows the occurrence rate of grid cells with a given temperature at a given height. An LTE equation of state results in three such horizontal bands, a non-equilibrium hydrogen ionization equation of state results in two, and a non-equilibrium hydrogen and helium ionization equation of state results in no horizontal bands. LTE ionization is described by a set of Saha equations. These equations give the ratio between two successive stages of ionization, and they are quite sensitive to temperature. According to these equations, a large change in the ionization state is associated with a minor change in the temperature. In the chromosphere, where the gas is partially ionized, the LTE ionization state will adjust to heating or cooling taking place, while the temperature remains roughly constant. The dark bands of figure 4.1 correspond to these "constant" temperatures. Further details on the implementation of non-equilibrium
Chapter 4. PhD project summary

Figure 4.1: Occurrence rate of grid cells with a given temperature and at a specific height. The panels show the results from four different simulation runs. They differ in how the equation of state is modelled and this is indicated by the label. LTE: all elements are in LTE. HION and LYA-HION are both solving the time dependent rate equations for hydrogen but with and without the Lyman-α transition. HELIUM: the time dependent rate equations for hydrogen and helium are solved. The simulation run with the LTE equation of state features three horizontal plateaus. The plateaus occur at the temperatures where ionization occur in LTE. The plateau corresponding to the ionization of hydrogen disappear in the simulation runs with a non-equilibrium hydrogen equation of state. The plateau featuring non-equilibrium of hydrogen and helium does not have plateaus.

helium ionization and its effects on the solar atmosphere can be found in Golding et al. (2016) which is given in chapter 6.

4.3 Formation of helium lines

Explaining spatial structure of 10830

The opacity of He I λ10830 in the chromosphere is known to be sensitive to the amount of impinging EUV photons emitted from the transition region and corona (see section 3.2). Solar images taken in the line show small scale structure. This suggests that localized EUV sources are at play, contributing to the formation of such structures. We investigate this by analysing a snapshot from a Bifrost simulation where the effects of non-equilibrium ionization of both hydrogen and helium are taken into account. The analysis is based on the radiative transfer calculations assuming statistical equilib-
4.3. Formation of helium lines

The formation of helium lines is carried out with the radiative transfer code Multi3d (Leenaarts & Carlsson, 2009). The emissivity due to EUV lines of the transition region and corona is self-consistently included in the calculations by assuming ionization equilibrium for the EUV emitting ions.

The analysis shows that the $\lambda 10830$ line deepens in regions that are experiencing strong EUV irradiation. In particular, chromospheric structures that are surrounded by a hot and dense transition region will show up as structures in the images. Further details about the study are found in Leenaarts et al. (2016) which is given in chapter 7.

On the formation of the EUV line intensity

Which mechanisms are dominating the formation of helium EUV resonance lines have been debated over the last decades (see section 3.1). To investigate possible mechanisms we use the new functionality of Bifrost to simulate the solar atmosphere including the effects of non-equilibrium ionization of hydrogen and helium. With a snapshot from the simulation we compute the emergent intensity in the helium EUV transitions and investigate the line formation in the model. The radiation transfer calculations are done with a version of Multi3d (Leenaarts & Carlsson, 2009) where we use the non-equilibrium helium ionization fractions from the model snapshot as constraints. With this code we obtain a solution of the radiative transfer equation that includes effects non-equilibrium ionization. The investigation shows that the enhanced intensity of He I $\lambda 584$ can be attributed to thick radiative transfer, and the enhanced intensity of the He II $\lambda 304$ and $\lambda 256$ lines can be attributed to non-equilibrium ionization and thick radiative transfer. Further details about the study are given in chapter 8.
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